# The origin of stars

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# Chapter 1

# Introduction

When we look up into the sky at night, the first thing we notice is a beautiful blanket of stars between all the blackness of space. Gazing into infinity, one of the first questions that must come to mind is where these bright dots that we call stars come from. This has to be the single most important reason that has propelled man to do astronomy. After millenia of research in physics and astronomy, one would expect that we have unraveled all the mysteries of this majestic puzzle. However, we have not. Like the vastness of the Universe, understanding the formation of stars is a complex beast by it self. We have made some progress over the years, of course, and our knowledge is gaining rapidly, still, we are like a child trying to understand the workings of a complicated toy. The study of star formation is therefore one of the longest standing research fields in astronomy, and dare I say, one of the most important as well.

Newly born stars form out of gas and dust from the remnants of their ancestors that enriched their environments during their lifetimes. Figure 1.1 shows one of the most famous star forming region images of modern times. Most of the star matter that is ejected into the space between the stars, the ISM (interstellar medium), comes from supernova explosions. Only massive stars, with masses above 8 solar masses, have the prerogative to perform this task. The majority of the heavy elements can only be formed and distributed in this way. This while the stars that have masses similar to, or lower than, our Sun are doomed to live for billions of years. Massive stars are therefore the heavy element factories of our Universe. Because of the continued evolutionary cycle, the initial conditions for each generation of stars will be different. Intuitively, this must have a significant influence on the early stages of star formation.

It is important to know how massive a star becomes when it forms. Or, what fraction of stars in a star forming region is massive enough to have a significant impact on its environment. This because they play key roles in many astronomical fields. Massive stars inject energy into the ISM through radiation, outflows, winds, and supernovae (Hensler 1999; Chappell & Scalo 2001). Knowing this, for example, will help us to better understand the star formation history of galaxies, the chemical enrichment of the early Universe, or even to predict the brown dwarf population in a cluster, and much more. It is therefore of great importance to quantify the relative numbers of stars in different mass ranges. This distribution of stellar masses at the moment of their creation is called the stellar initial mass function.

This thesis tries to give insight into the evolution of interstellar clouds and its dependence on the initial and ambient environmental conditions in the most extreme environments of our Universe. It looks into the fragmentation properties of molecular clouds and into the formation of stars using numerical simulations, thereby focussing on its dependence on metallicity, rotation, and feedback effects (mechanical and radiative). Trying to answer the universality conundrum of the initial mass function in extreme environments is a key aspect and the main goal of this work.

**Figure 1.1:** Pillars of Creation. The giant pillars consist of molecular gas and dust, which are so dense that the interior gas contracts gravitationally to form stars. Intense UV radiation from nearby massive stars evaporates the surfaces of these clouds to give them their present shape. This image, which lies inside the Eagle nebula at a distance of 2 kpc away from us, was taken with the Hubble space telescope in 1995. Image credit: NASA, ESA, STScI, HST, J. Hester and P. Scowen (Arizona State University).

## 1.1 The theory of star formation

Star formation is a complex dynamical process. Stars are known to form inside huge gas clouds that are rich in various kinds of molecules. It starts out with clumps of gas and dust (if present) merging with each other and held together by their small gravitational forces. When enough mass piles up, the gravitational pull increases and the cloud starts to contract. There is, however, pressure which prevents a cloud to collapse under its own gravity. Not only thermal and radiation pressure, but also magnetic fields and bulk motions are sources that can prevent collapse. Only when gravity dominates all these counter forces, collapse will occur and continue to form a star. An image to this effect is shown in Fig. 1.2.

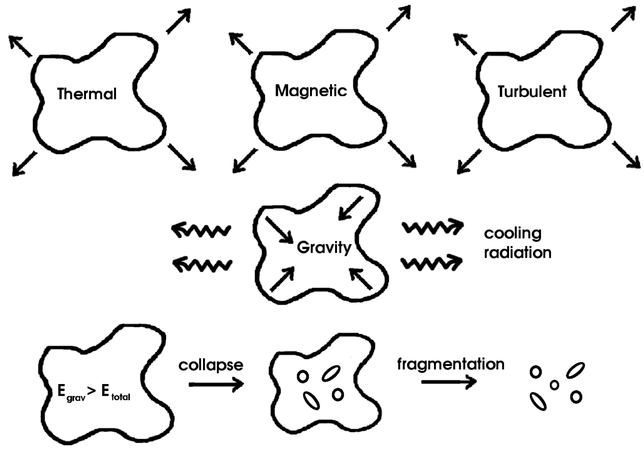


**Figure 1.2:** The collapse of a molecular cloud. Gravitational pull needs to be stronger than all the other forces in order to initiate collapse and subsequent fragmentation.

One can calculate how much (minimal) mass is needed for gravity to dominate all other forces and start the formation process by using the Jeans mass formula

$$M_J = \left(\frac{3}{4\pi}\right)^{\frac{1}{2}} \left(\frac{5k_B}{G\mu m_H}\right)^{\frac{3}{2}} \frac{T^{\frac{3}{2}}}{\rho^{\frac{1}{2}}}, \tag{1.1}$$

where, $\mu$ is the mean molecular weight, $m_H$ is the mass of a hydrogen atom, G is the gravitational constant, $k_B$ is the Boltzmann constant, T is the gas temperature, and $\rho$ is the mass density. As can be seen from this equation, the critical mass is only a function of temperature and density. Increasing density and decreasing temperature lead to lower Jeans masses which, in turn, will lead to fragmentation into smaller cores. It is possible to formulate this equation in a more practical form as given by Frieswijk et al. (2007), that is,

$$M_J \simeq \frac{90}{\mu^2} M_\odot \left(\frac{n}{1\,\mathrm{cm}^{-3}}\right)^{-\frac{1}{2}} \left(\frac{T}{1\,\mathrm{K}}\right)^{\frac{3}{2}}. \tag{1.2}$$

When an interstellar cloud of gas is in hydrostatic equilibrium, the internal energy is in balance with the potential energy of the self-gravitational force. The virial theorem states that twice the average total kinetic energy must equal n times the average total potential energy for a system to be in a stable statistical equilibrium. If the system consists of matter held together by its own gravity, then, n equals -1. From the virial theorem, one can simply derive the Jeans equation. If we assume that the kinetic energy, $E_k$, is dominated by the thermal energy, and the potential energy, $E_p$, by the gravitational energy, then the virial theorem

$$2E_k = -E_p \tag{1.3}$$

can be formulated as

$$f_1 N k_B T = f_2 \frac{GM^2}{r}, \tag{1.4}$$

where $f_1$ is a constant that depends on the degrees of freedom of the system, which is 3/2 for a monatomic gas, and $f_2$ is a constant which depends on the shape of the body and is 5/3 for a sphere. N is the total number of particles, which can be substituted by $M/\mu m_H$ in Eq. 1.4 and r, the radius of the cloud, can be substituted by $(3M/4\pi\rho)^{1/3}$, such that Eq. 1.1 is recovered.

It must be noted that, in these equations, the assumption is made that the system constitutes an ideal, isothermal gas with negligible magnetic field effects and negligible bulk motions. While r is taken as the average radius of the cloud and where the density does not depend on the radius. We can eliminate some of these assumptions and can obtain the Jeans mass again by expressing the Jeans length, $\lambda_J$, in terms of sound speed $c_s$, where $c_s^2 = \gamma P/\rho$ and $c_s^2 = \gamma T$ in an ideal gas. Here, P denotes the pressure and $\gamma$ is a factor that depends on the equation of state, which will be elaborated on in section 1.3. Using these expressions, we can now start to derive the magnetic and the turbulent Jeans masses. The Jeans length is given by

$$\lambda_J = \left(\frac{\pi c_s^2}{G\rho}\right)^{\frac{1}{2}}. \tag{1.5}$$

For a pressure of the form: $P = k_B \rho T/\mu m_H$, Eq. 1.1 is reobtained. However, if magnetic fields were to dominate the pressure, where $P_{mag} = B^2/8\pi$ and thus $c_s^2 = \gamma B^2/8\pi\rho$, the Jeans mass would become

$$\begin{aligned} M_J^B &\propto \lambda_J^3 \rho = \left(\frac{\pi c_{s,B}^2}{G}\right)^{\frac{3}{2}} \rho^{-\frac{1}{2}} \\ &\simeq \frac{1.24 \times 10^7}{\mu^2 \gamma^{-\frac{3}{2}}} M_\odot \left(\frac{n}{1\ \mathrm{cm}^{-3}}\right)^{-2} \left(\frac{B}{3\mu G}\right)^3, \end{aligned} \tag{1.6}$$

where $B = 3\mu G^*$ is the typical magnetic strength in the ISM (Subramanian & Barrow 1998; Schleicher et al. 2009).

The observed turbulent motions in cloud cores certainly contribute as a counterforce to the gravitational pull. If the turbulence is isotropic, then the Jeans mass can simply be computed using an effective sound speed $c_{s,eff}^2 = c_s^2 + \frac{1}{3} v_{rms}^2$. In the case that turbulence dominates, the Jeans mass then becomes (Larson 1998; Elmegreen 1999; Spaans & Silk 2005; Frieswijk 2008)

*Here, $\mu G$ stands for micro Gauss. Not to confuse with grav. constant G and mean molecular mass $\mu$.

$$
\begin{aligned}
M_J^{\rm turb} &\propto \lambda_J^3 \rho = \left(\frac{\pi c_{\rm s,turb}^2}{G}\right)^{\frac{3}{2}} \rho^{-\frac{1}{2}} \\
&\simeq \frac{2.41 \times 10^4}{\mu^{\frac{1}{2}}} M_\odot \left(\frac{n}{1\ {\rm cm}^{-3}}\right)^{-\frac{1}{2}} \left(\frac{v_{\rm rms}}{1\ {\rm km/s}}\right)^3 . \qquad (1.7)
\end{aligned}
$$

When each of these conditions is fulfilled and collapse has started, one or more stars can be formed. It is, of course, not as simple as that. There are many other processes that the system goes through before reaching the final stage of a star. During collapse, angular momentum will increase, the density and pressure will rise, shocks will occur, energies will interchange, internal heat can be radiated away or trapped, and the molecular cloud might fragment. Feedback effects, such as, chemical, radiative, or mechanical, from either internal or external sources, will also play key roles. It is in these processes that the stage is set for the eventual masses of stars.

## 1.2 The interstellar medium

The interstellar medium is the medium between the stars inside galaxies, and interstellar space is not empty. The ISM is filled with pockets of gas with relatively low density ($\sim 1$ cm$^{-3}$), on average, and is mixed with dust. By comparison, the air on earth has a number density of $2.5\times10^{19}$ cm$^{-3}$. This means that the average human lung can inhale more atoms per hour than the average amount of interstellar gas in a volume half the size of the moon. As such, one needs gigantic gas clouds to be able to make even a few stars. There are such clouds. Typical sizes for molecular clouds span a range from 0.1 pc (stellar cores) to 100 pc (giant molecular clouds, Frieswijk 2008).

The primary sites for star formation are molecular clouds. These are thought to form out of the remnants of supernova explosions, outflows and winds from stars, and the gas reservoirs in the interstellar and the intergalactic medium (Ballesteros-Paredes et al. 1999; Hartmann et al. 2001; Klessen et al. 2005a; Heitsch & Hartmann 2008; Dobbs & Bonnell 2008; Dobbs 2008). A molecular cloud comprises atoms, molecules, ions, and dust which make it a rich chemical system. Molecules form in the gas phase or on dust grains and allow low temperatures to be reached (Omukai et al. 2005; Cazaux et al. 2005a; Cazaux & Spaans 2009; Dulieu et al. 2009). The composition of these clouds depend on the ambient conditions like temperature, pressure, density, and metallicity. Multiple phases can exist within one cloud with different species and temperatures.

Molecular clouds in the Universe have different shapes and sizes. Fig. 1.3 depicts three very well known gaseous nebulae. They are morphologically quite different from one another. The density is also not constant throughout interstellar clouds. Hot ionized regions usually occupy the low density medium $n \ll 1$ cm$^{-3}$, while molecular cores reside in rather dense locations $n \geq 10^4$ cm$^{-3}$.

**Figure 1.3:** Three interstellar clouds. On the top, the Orion nebula is shown. On the bottom left is the Tarantula nebula and on the bottom right is the Rosette nebula. Image credit: ESA, Herschel, NASA.

The density profile is not known for the larger molecular clouds ($r \gtrsim 0.1$ pc). They have bulky shapes with a perturbed density structure. Besides, it is not likely that a well-developed density profile exists before virialization. Modelers working in this field commonly assume either a flat density profile or some random perturbation, while providing the models with realistic velocity profiles, which naturally evolve into the observed sub-structures. According to some astronomers, the distribution of the dense regions in such clouds could be the precursor of the stellar initial mass function (Motte et al. 1998; Motte & André 2001; Goodwin et al. 2008; Simpson et al. 2008; André et al. 2010), and perhaps the stellar IMF is directly linked to its dense core mass function, as it seems for the Pipe nebula (Lada et al. 2008). The dense cores ($r \sim 0.1$ pc), on the other hand, do have a well defined density structure. Observational evidence shows that pre-stellar dense cores increase their density towards the center and then reach a plateau. Commonly used density profiles when modelling these cores are either a $\rho \propto r^{-2}$ profile or a Plummer-like, $\rho = \rho_0 \left(1 + r^2/a^2\right)^{-\eta}$, profile, where $\eta = 5/2$ is the classical Plummer sphere and, a is a scale parameter which sets the size of the core.

### 1.2.1 Fragmentation

A large cloud that is collapsing has the tendency to fragment. Gas above the critical mass ($M > M_J$) will go into a free-fall when there is no significant counter pressure. This is likely, at least initially, since molecular clouds are optically thin so that radiation can easily escape, cooling the cloud and releasing any pressure. The free-fall time only depends on the gas density as follows

$$\tau_{ff} = \sqrt{3\pi/32G\rho}. \tag{1.8}$$

Each supercritical overdensity will have a different free-fall time. The overdensities will therefore contract at different rates. As the molecular cloud collapses, it will break into smaller and smaller pieces in a hierarchical manner, until the fragments reach their Jeans mass. As the density increases, each of these fragments will become increasingly more opaque and are thus less efficient at radiating away the gravitational potential energy. This raises the temperature of the cloud and inhibits further fragmentation.

The fragments might normally merge with one another if the increasing angular momentum did not prevent this. As the cloud shrinks due to gravity, it spins faster. This hinders the collapse and favors fragmentation of the cloud. The main mechanism that causes molecular clouds to have overdensities in the first place as well as sustaining an asymmetric velocity structure is thought to be turbulence. The interplay between gravity and turbulence which enhances fragmentation is also described as gravoturbulent fragmentation (Klessen & Ballesteros-Paredes 2004).

### 1.2.2 Turbulence

Molecular clouds are observed to be quite turbulent (Larson 1981; Falgarone et al. 2001; Caselli et al. 2002; Heyer & Brunt 2004; Brunt et al. 2009; Brandenburg & Nordlund 2011). It is not precisely known where the interstellar turbulence comes from, however, there are some reasonable ideas. Turbulence can originate from supernova explosions which send shock waves into the ISM at very high speeds. When these shock waves come in contact with an object, or another shock wave, they can dissipate and cause turbulent motions. This is thought to be one of the main drivers of turbulence. Another culprit that can incite turbulence is infalling matter from the intergalactic medium (IGM). Gaseous matter can flow into the ISM from the IGM due to the potential wells of galaxies. Upon mixing with the interstellar matter, large scale turbulence is created. Recent discoveries show that in order to maintain long term star formation, continues gas supply is needed. Huge reservoirs of gas are available in the IGM, which is the most logical source of this fuel (Bauermeister et al. 2010). Other astrophysical processes acting on large scales, including magneto-rotational instability, or spiral shock forcing, are also thought to be viable candidates for the generation and maintenance of molecular cloud turbulence (Brunt et al. 2009). Gravitational instabilities during contraction are strong drivers of turbulence as well, however, they act on smaller scales and could not instigate the initial turbulence. Fig. 1.4 shows a picture of a simulated turbulent cloud.

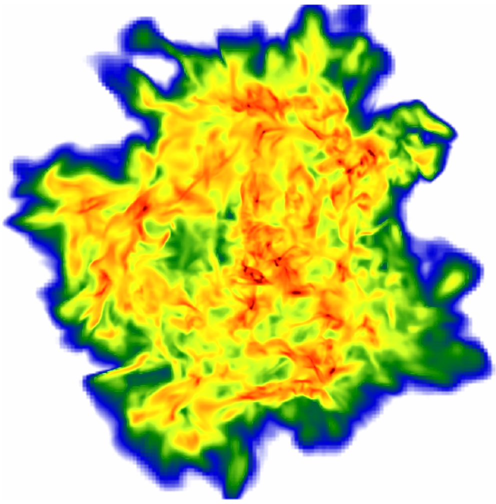

**Figure 1.4:** A simulated turbulent molecular cloud. Image by Hocuk & Spaans (2010a).

The velocity scaling of turbulent motions appears to be similar in regions with varying intensity of star formation (Brunt & Heyer 2002a,b). This indicates that the velocity scaling is inertial, and driven mostly by energy input at large scales, rather than local driving by on-going star formation (Brandenburg & Nordlund 2011). To this end, it is thought that turbulence caused by stellar feedback, e.g., radiative or mechanical, although an integral part of the formation processes, is not strong enough to drive the large scale turbulence in molecular clouds (Brunt et al. 2009). This while jets from stellar outflows might produce strong shocks or disturbances which can trigger or disrupt local gravitational collapse. All of these mechanisms do contribute a lot of mechanical energy into the interstellar medium and intensify the existing turbulence.

The turbulence in interstellar clouds is found to behave like a power-law (Larson 1979, 1981; Ossenkopf & Mac Low 2002), where the largest scales contain the highest energies. The velocity scaling is of the form

$$(\sigma_{\rm v} =) \triangle {\rm v} \propto \ell^m, \tag{1.9}$$

where $\sigma_{\rm v}$ and $\triangle$v denote the velocity dispersion, $\ell$ the length scale, and $m$ the scaling power. The value of $m$ for compressible fluids would be $m = 1/2$ in this notation, while incompressible, Kolmogorov turbulence has a value of $m = 1/3$. The power spectrum commonly used in numerical simulations directly follows from this velocity scaling. The power spectrum of turbulent motions as observed in molecular clouds (Larson 1981; Myers & Gammie 1999; Heyer & Brunt 2004) scales as

$$\rm P(k) \propto k^{-4}, \tag{1.10}$$

where $\rm k = 2\pi/\ell$ is the wavenumber.

It is sometimes preferred to write the power spectrum in the form of an energy relation, also called the energy spectrum. Considering that the relation between the energy spectrum and the power spectrum is $E(k) = 2\pi k^2 P(k)$, the energy spectrum scales with the wavenumber as $E(k) \propto k^{-2}$.

### 1.2.3 Metallicity and Dust

Metallicity (Z) is a term used by astronomers and describes the relative abundances of elements. All elements heavier than helium are collectively labeled as metals. Metallicity increases as the ratio of heavy elements (metals) with respect to helium and hydrogen becomes larger. The metallicity of interstellar clouds in the solar neighbourhood is similar to that of our Sun ($Z_\odot$). Of the gas in the ISM, 93% of the atoms (by number) are hydrogen and 7% are helium, with $< 1\%$ of atoms being elements heavier than hydrogen or helium. Metallicity is an important factor in star formation, since the creation and propagation of radiation through a medium is affected by absorption, emission, and scattering processes, which, in turn, are all dependent on the composition of the medium.

All species (atoms, molecules, and ions) can absorb or emit radiation through various means (free-free, bound-free, or bound-bound processes). Generally, species with a higher atomic number, and with more degrees of freedom, have a wider range of possible transitional states. Therefore, they are more efficient in releasing the internal energy of a system. A high metallicity content of a molecular cloud will allow it to cool much faster and allow lower temperatures to be reached. Until local thermodynamic equilibrium (LTE) is obtained, the rate of cooling scales with the square of the density. The cooling rate depends on a cooling function as follows

$$\Lambda_{cr} = 4\Lambda_{cf} n_p n_e \underbrace{\left(\frac{n_c}{n_c + n_e}\right)}_{\text{correction for LTE effects}}, \tag{1.11}$$

where $n_p$ is the number density of protons, $n_e$ is the number density of electrons, $n_c$ is the critical density around which the transition to LTE occurs, which is $\sim 10^3$ cm$^{-3}$, and $\Lambda_{cf}$ is the cooling function. A cooling function depends on various quantities, e.g., on the atomic and molecular composition of the gas, on the properties of dust, and on physical conditions like temperature and density of the medium. Fig. 1.5 shows four cooling functions for different metallicities created with the Meijerink & Spaans (2005) code.

Cooling is only important in the evolution of the molecular cloud if thermodynamic processes occur on a timescale shorter than the collapse timescale (Eq 1.8). The cooling timescale is given by

$$\tau_{cool} = \frac{3k_B \rho T}{2m_H \Lambda_{cr}}. \tag{1.12}$$

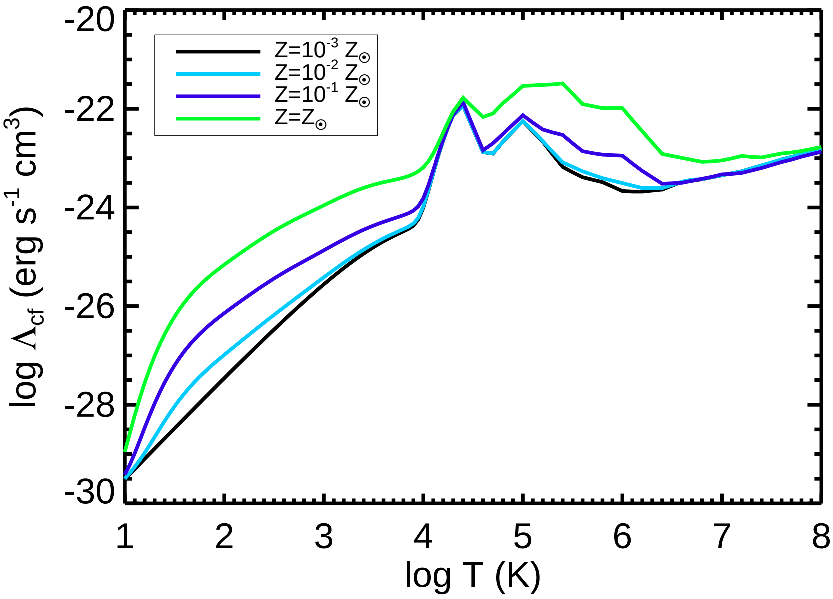


**Figure 1.5:** Cooling functions. Four detailed functions created by Meijerink & Spaans (2005) corresponding to four metallicities $Z = Z_{\odot}$, $10^{-1}$ $Z_{\odot}$, $10^{-2}$ $Z_{\odot}$, and $10^{-3}$ $Z_{\odot}$.

Aside from alleviating a molecular cloud from its internal energy by radiating it away, metals absorb radiation much easier as well. They cause a cloud to reach higher opacities much quicker during collapse. This will result in the trapping of radiation and the molecular cloud will no longer be able to cool down further, thereby increasing its temperature and pressure as it contracts. Gas becomes optically thick at number densities of $\sim 10^{16}$ $cm^{-3}$, if it is fully atomic, while it becomes optically thick around $10^{12}$ $cm^{-3}$ when it has solar metallicity (Machida et al. 2009). For a collapsing molecular cloud, the interplay between heating and cooling is vital in determining the state at which stars form (Hocuk & Spaans 2010b).

Dust also plays an important role in the formation of molecules that affect the star formation processes (Cazaux et al. 2010, 2011a,b). Molecular clouds contain about 1% dust, by mass (Cazaux 2004). In astronomy, dust is a term used for aggregates of material, like chains of molecules, such as PAHs (poly-aromatic hydrocarbons) and small grains, held together by various bonds. In order to form dust grains one needs atoms with unique bonding properties, like carbon or silicon, because it is easy to form chains with them. It is therefore obvious that the dust content must depend on the metallicity of the medium and is thought to be tightly correlated with it (Cazaux & Spaans 2009). In Fig. 1.6, an image of such a grain is shown.

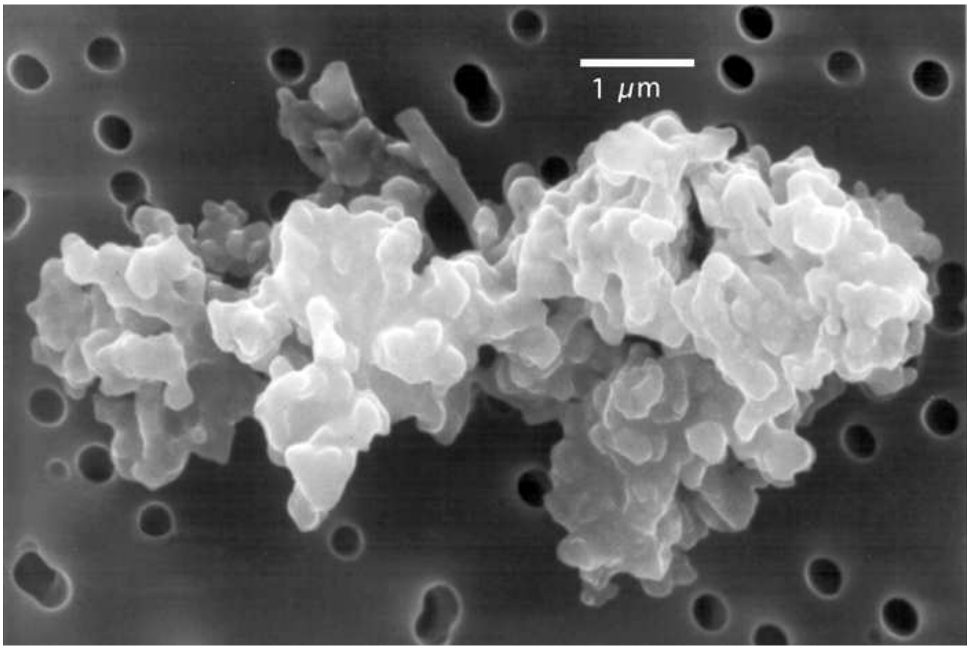


**Figure 1.6:** A porous chondrite dust particle. Courtesy of E.K. Jessberger.

Dust will also affect the thermal balance of the gas. When gas and dust are collisionally coupled ($n_H \gtrsim 10^{4.5}$ cm$^{-3}$ for solar metallicity), gas can either cool down or heat up on the dust grains. More importantly, dust can enhance the formation of molecules, especially $H_2$, by acting as a catalyst at temperatures above $T > 20$ K (Cazaux et al. 2005b). At lower temperatures, gas can freeze onto the dust grains, thereby locking the species up in them until it is warm enough again for evaporation. This regulates the temperature of a system until a stable point is reached. Despite the fact that the metallicity of the system remains unchanged, creating new and more complicated species helps to cool the system much more efficiently.

## 1.3 Equation of state

The equation of state (EOS) describes the relation of the state variables in a system to one another. In fluid dynamics, typically, the relation of pressure against density and temperature is considered. The pressure of an ideal gas scales linearly with temperature and density. More precisely, it behaves in the following manner

$$P = \frac{N_A k_B \rho T}{\mu}, \tag{1.13}$$

where $N_A$ is Avogadro's constant, $k_B$ is the Boltzmann constant, T is the gas temperature, and $\mu$ is the average mass per particle.

It is also possible to relate pressure and density independently from temperature. This is by assuming a polytropic equation of state. A polytropic process is a thermodynamic process that obeys the relation

$$PV^n = \text{Constant}, \tag{1.14}$$

where n is the polytropic index and V is the volume. But, how do we get to this equation? Sometimes, it is possible to write down the equations which describe a physical system, but for which solutions cannot be derived analytically. In many cases, the only way to solve a specific problem will be using numerical, computer based, methods. In the case of a self-gravitating, spherically symmetric fluid, however, it is possible to analytically solve the equation characterizing the system, also known as the Lane-Emden equation, by using a polytrope. For example, if we take the equation for hydrostatic support in terms of the radius variable r

$$\frac{dP(r)}{dr} = -\frac{GM(r)\rho(r)}{r^2}, \tag{1.15}$$

in which G is the gravitational constant and M(r) is the mass at radius r, and take the derivative with respect to r, we get

$$\frac{d}{dr}\left(\frac{r^2}{\rho}\frac{dP}{dr}\right) = -G\frac{dM(r)}{dr}. \tag{1.16}$$

Now, if we solve the right hand side, we obtain

$$\frac{1}{r^2}\frac{d}{dr}\left(\frac{r^2}{\rho}\frac{dP}{dr}\right) = -4\pi G\rho. \tag{1.17}$$

This second order differential equation can be solved using a polytropic EOS of the form

$$P \propto \rho^{1+1/n} = \rho^{\gamma}, \tag{1.18}$$

with $\gamma$ being the polytropic exponent.

A polytropic EOS is not really independent of temperature. The temperature dependency is rather hidden in the exponent $\gamma$, which is an important parameter that explains the behaviour of the system. $\gamma$ is typically considered to be a function of physical parameters or ambient conditions, such as, gas temperature, dust temperature, radiation intensity, velocity field, chemical composition, metallicity, and magnetic field (Spaans & Silk 2000). A $\gamma = 1$ would, for example, tell us that the system is isothermal, since the EOS becomes $P \propto \rho$ now. Whereas $\gamma = 5/3$ describes an adiabatic system, typical for high density gas and inside main sequence stars. When we consider Eqs. 1.13 and 1.18, take their log and then the derivative, i.e.,

$$\begin{aligned} \mathrm{dlogP} &= \mathrm{dlog}\rho + \mathrm{dlogT} \\ \mathrm{dlogP} &= \gamma \mathrm{dlog}\rho, \end{aligned} \tag{1.19}$$

we acquire the relation for $\gamma$,

$$\gamma = 1 + \mathrm{dlogT}/\mathrm{dlog}\rho. \tag{1.20}$$

This tells us that $\gamma$ depends on the details of heating and cooling of a system through these derivatives and, as such, depends implicitly on the radiative transfer effects and the changes in the chemical composition. This description of the EOS is true as long as the heating and cooling terms in the energy equation balance out on a shorter timescale than the gas dynamics do (Scalo & Biswas 2002; Spaans & Silk 2005).

One interesting feature of the polytropic EOS is the $\gamma$ at which the Jeans mass remains constant. Since the Jeans mass, $M_J$, scales as $T^{\frac{3}{2}}/\rho^{\frac{1}{2}}$, and one can rewrite the temperature as a function of density by combining Eqs. 1.13 and 1.18, that is, $T = \rho^{\gamma-1}$, the Jeans mass then becomes proportional to

$$M_J \propto \rho^{\frac{3}{2}\gamma-2}. \tag{1.21}$$

This means that the Jeans mass becomes independent of the variables and remains constant when $3/2\gamma - 2 = 0$. A $\gamma$ of 4/3 fulfills this condition. One can see from this that for a $\gamma$ lower than 1.33, the Jeans mass decreases as the gas contracts and the density increases. Thermal pressure can no longer stop the collapse and a molecular cloud in such a state will be highly susceptible to fragmentation.

A value of $\gamma$ below 1, but still greater than 0, would mean that the temperature decreases as the density increases, while the pressure still increases at a slow rate. This can only occur in a system that is losing its internal energy by any means, like radiating it away. This is frequently seen in optically thin interstellar clouds. As such, the softness of $\gamma$ plays a major role, at a very early stage, in the fragmentation properties and the evolution of molecular clouds (Spaans & Silk 2000; Li et al. 2003; Klessen et al. 2005b; Jappsen et al. 2005). A negative $\gamma$ is an unlikely state in which even the pressure decreases with increasing density. Such a state cannot be sustained for long periods of time, but is known to occur in explosions.

## 1.4 The initial mass function

The initial mass function (IMF) is a distribution of stellar masses versus their number in a given volume of space. It is an empirical function that is observed to behave like a power-law for masses above a few tenths of a solar mass. This renowned function is obtained through a logarithmically binned histogram of the initial masses of stars. Fig. 1.7 shows an example plot of an IMF that is representative of the star forming conditions in our solar neighbourhood.

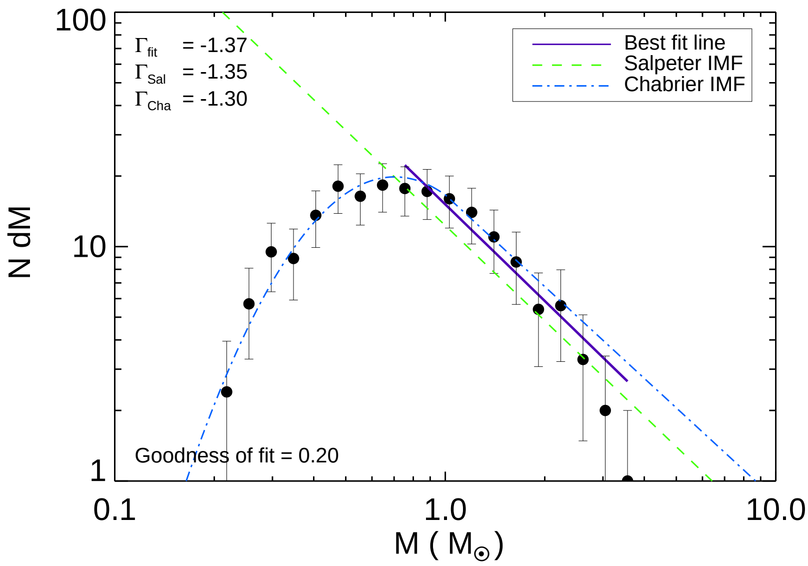


**Figure 1.7:** A simulated IMF of the Milky Way. This mass function is obtained from numerical simulation using conditions of the ISM similar to our Galaxy. Two commonly used IMFs, the Salpeter IMF and the Chabrier IMF, are overplotted in this figure to highlight the good match.

The IMF was first proposed by Edwin Salpeter in his famous paper (Salpeter 1955). He found from his observations of stars that plotting the number of stars against their mass resulted in a power-law. The functional form that he proposed for the IMF was the following

$$\frac{dN}{dM} \propto M^{-\alpha}, \tag{1.22}$$

with N the number of stars in a mass range dM, $\alpha$ the power-law index above the characteristic mass of $\sim 0.3\ M_\odot$, also known as the turn-over mass. Edwin Salpeter empirically found that the power-law index has a value of $\alpha = 2.35$. At that time, observational constraints made it difficult to resolve stellar masses below the mass of our Sun and because of this, he could not see a turn-down of the mass function at lower masses. Until today, the value of the power-law index has remained remarkably unchanged above $1\ M_\odot$ and is now known as the Salpeter slope.

It is often more practical to formulate the IMF in a logarithmic form. The following steps show the derivation to this end. To get the number of stars within the mass range M+dM, one can integrate Eq. 1.22 as follows

$$N = \int dN \propto \int M^{-\alpha}\, dM = (-\alpha+1)^{-1}\, M^{-\alpha+1}. \tag{1.23}$$

Now, by taking the logarithm on each side and looking only at the change as a function of (log) mass, thereby losing the constants, one will end up with the following equation;

$$\frac{\mathrm{dlogN}}{\mathrm{dlogM}} = -\alpha + 1 = \Gamma. \tag{1.24}$$

Here, $-\alpha + 1$ is replaced by $\Gamma$, in which $\Gamma$ defines the slope above the characteristic mass. This makes the slope of the Salpeter function $\Gamma = -1.35$. Eq. 1.24 is the generally used convention by many astronomers and the adopted form throughout this thesis.

In the following figure (Fig. 1.8), we illustrate three well known IMFs. These are, the Salpeter (1955) IMF (purple), the Chabrier (2003) IMF (red), and the Kroupa (2001) IMF (green). While the Salpeter IMF has a slope of $\Gamma = -1.35$, the Chabrier IMF is a log-normal function up to 1 solar mass. Above 1 solar mass, it follows a power-law with a slope of $\Gamma = -1.3$. The Kroupa IMF on the other hand, is a canonical IMF with three different power-law slopes. These are: $\Gamma = 0.7$ between 0.01 $M_\odot$ and 0.08 $M_\odot$, $\Gamma = -0.3$ between 0.08 $M_\odot$ and 0.5 $M_\odot$, and $\Gamma = 1.3$ between 0.5 $M_\odot$ and 50 $M_\odot$.

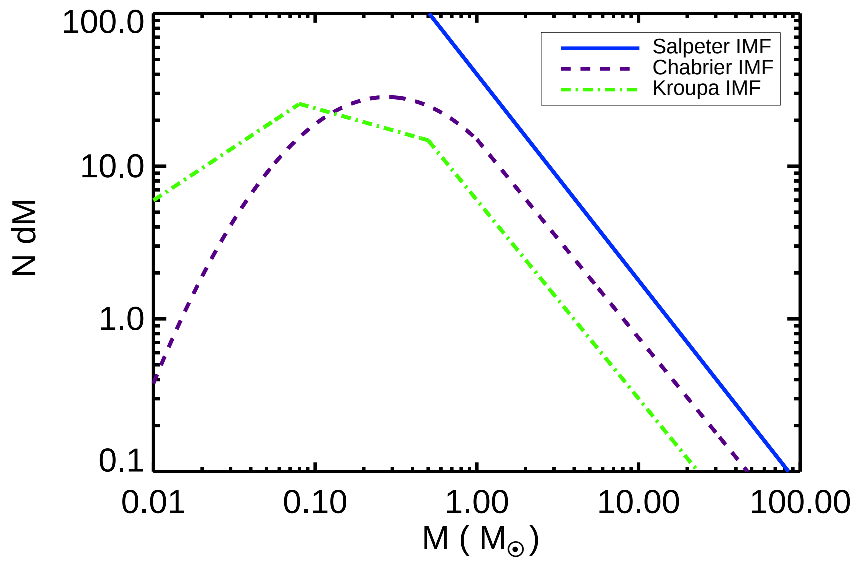


**Figure 1.8:** Three different initial mass functions. The purple line shows the Salpeter IMF. The red dashed line shows the Chabrier IMF. The green dot-dashed line shows the Kroupa IMF. The Y-axis is in arbitrary units. Image created by author.

The initial mass function offers the great benefit of predicting the likelyhood of stellar masses at the time of their formation. It has become an important diagnostic tool and is of fundamental importance for many astronomical fields. The IMF is used in areas, such as, galaxy formation and evolution, studies of chemical enrichment, the prediction of brown dwarfs, energetic feedback into the ISM, and in just about anything that studies large quantities of stars (Zoccali et al. 2000; Kroupa 2002; Bonnell et al. 2007). It is therefore appropriate to state that one of the main goals for a theory of star formation is to understand the origin of the stellar initial mass function.

It has been theorized that this distribution is universal in nature, nicely following a Salpeter slope (Salpeter 1955). Numerous observations of the solar neighbourhood support this hypothesis (Chabrier 2003; Sabbi et al. 2007; Elmegreen et al. 2008) and has gained a lot of interest since this idea first arose. The shape of the distribution has been further refined by other astronomers, especially improving upon the lower mass end of the function (Miller & Scalo 1979; Kroupa 2001; Chabrier 2003). A universal IMF would be quite useful, as one can imagine, since it would eliminate one major uncertainty in a lot of astronomical fields.

As charming as the idea is, the story does not end there. Several recent studies, of mainly extragalactic objects, have started to show us the other side of the story. At the turn of the century, astronomers were beginning to notice deviations in their mass functions from a standard Salpeter shape (Baugh et al. 2005; Nayakshin et al. 2007; Parra et al. 2007; Davé 2008; Wilkins et al. 2008; van Dokkum 2008; Elmegreen 2009). These came from measurements of abundance patterns in extragalactic bulges (Ballero et al. 2007, 2008), enhancement of far infra-red luminosities in interacting galaxy systems (Brassington et al. 2007), mass-to-light ratios of ultra-compact dwarf galaxies (Dabringhausen et al. 2009a), NaI and FeH band spectra in luminous elliptical galaxies (van Dokkum & Conroy 2010, 2011), and many others. Our Galaxy center also shows signs of variations in the IMF (Figer 2005a,b; Paumard et al. 2006; Espinoza et al. 2009; Elmegreen 2009; Bartko et al. 2010). Complementary to these observations, many numerical studies support a non-universal IMF as well (Klessen et al. 2007; Hsu et al. 2010; Krumholz et al. 2010; Girichidis et al. 2011). Yet, they were much debated and solid evidence is still lacking. These developments in the last decade have driven astronomers towards a crossroads, and has divided them, which was the beginning of the IMF's universality conundrum.

Do all star forming events give rise to the same distribution of stellar masses? Is star formation essentially a self-regulating process or is fragmentation the process by which stellar masses are fixed? Certainly, the details of either process depend on the physical conditions of the cloud of gas and dust from which the stars form. The unanswered question is: how sensitively does the distribution of stellar masses depend on the initial conditions in the natal environment? Our ultimate goal is to understand the physical mechanisms that are responsible for the origin of stellar masses.

## 1.5 Extreme environments

The process of star formation is poorly understood in extreme environments. It is uncertain if stars form in the same way everywhere and if the IMF is similar to our Galaxy or that star formation is severely affected by the harsh, non-Milky Way ambient conditions. Our Universe harbors many regions which are quite extreme. An extreme environment in space exhibits conditions that are challenging to the formation of stars. These may include high temperatures, high pressures, strong radiation fields, powerful turbulence, strong gravity, or very low metallicities. Studying the evolution of gaseous clouds and star formation in extreme environments will help us understand the processes of star formation by determining its limits.

### 1.5.1 Active galactic nuclei

Active galaxies harbor supermassive black holes with masses of $\gtrsim 10^6$ $M_\odot$ in their centers. Active galactic nuclei (AGN) are the most luminous sources of electromagnetic radiation in the Universe. These accrete matter from the inner parsecs of galactic centers. Accretion rates can go up to their Eddington limit, if magnetic fields are not significant, but are typically $\sim$10% of this limit (Meijerink et al. 2007). It is also possible to surpass this limit in certain circumstances which might play an important role in the gas accretion process in host galaxies (Kawakatu & Ohsuga 2011). The Eddington rate is obtained when the hydrostatic equilibrium equation, as given in Eq. 1.15, is set equal to the continuum outward radiation pressure

$$\frac{dP_{rad}}{dr} = -\frac{\kappa\rho}{c}F_{rad} = -\frac{\sigma_T\rho}{m_p c}\frac{L}{4\pi r^2}, \tag{1.25}$$

where $\sigma_T = 6.65 \times 10^{-25}$ cm$^2$ is the Thomson scattering cross section for the electron and the gas is assumed to be purely made of ionized hydrogen, $\kappa$ is the opacity coefficient of the stellar material, $F_{rad}$ is the radiation flux, L is the luminosity, c is light speed, and $m_p$ is the proton mass. The Eddington luminosity then becomes

$$L_{Edd} = \frac{4\pi G M m_p c}{\sigma_T} \text{ erg s}^{-1}, \tag{1.26}$$

while the accretion rate can be written as

$$\dot{M} = \frac{dM}{dt} = \frac{2L_{Edd}}{c^2} = \frac{8\pi G M m_p}{c\sigma_T} \text{ g s}^{-1}. \tag{1.27}$$

The matter around black holes is heated up through the accretion process to millions of degrees. This heating leads to the emission of highly energetic radiation, such as UV photons (E > 5 eV) and X-rays (E > 1 keV). The accretion disks shine so brightly that they are visible from great distances, which renders them easy to spot with telescopes. Besides strong radiation fields, supermassive black holes impart a strong gravitational field as well. Molecular clouds in the inner 10 pc of active galaxies will be strongly affected by this gravitational pull as well as the irradiation by X-rays and UV (Meijerink et al. 2007; Hocuk & Spaans 2010a).

### 1.5.2 Starbursts

Starbursts are regions of space with an unusually high rate of star formation. Star formation rates (SFR) are observed as high as several hundreds to thousand solar masses per year and on the scales of a Galaxy (Sanders & Mirabel 1996; Smail et al. 1997; Hughes et al. 1998; Genzel et al. 1998). Massive stars are thought to commonly form in these places and an exceptional amount of UV radiation (E = $5 - 100$ eV) can be seen. The output of UV radiation is dominated by the O and B stars as only the hottest stars produce them. The supernova rate is also generally high in starbursting galaxies. This will in turn lead to an enhanced cosmic ray rate, since cosmic rays are thought to be produced mainly in supernova remnants (Papadopoulos et al. 2011). Figure 1.9 shows the famous starburst 'Antennae Galaxies'.

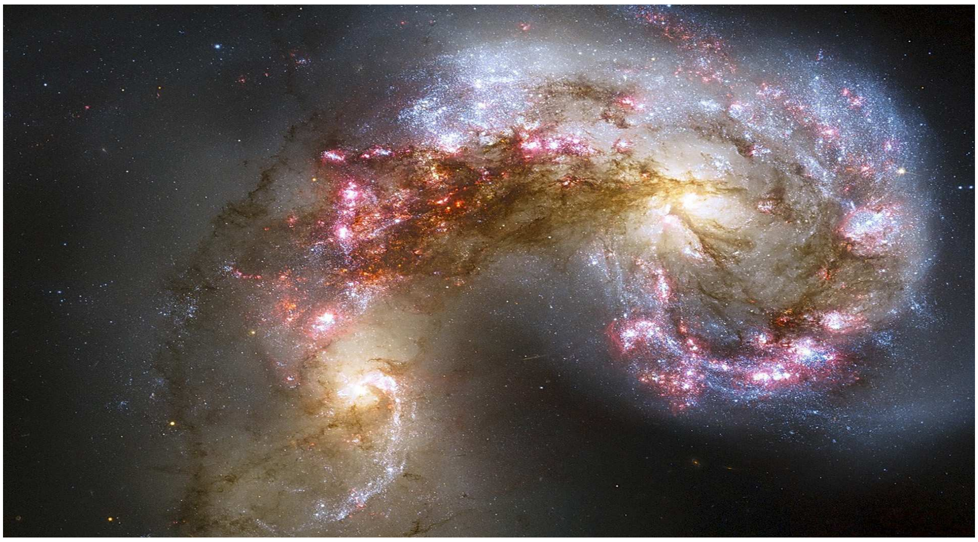

**Figure 1.9:** The Antennae Galaxies are an example of a very high starburst galaxy occurring from the collision of NGC 4038/NGC 4039. Credit: NASA/ESA

Over the years, mounting empirical evidence has been found that there is a correlation between nuclear activity and star formation over a wide range of redshifts (Kennicutt 1998; Trichas et al. 2009; Wild et al. 2010). It is thought that there is a tight relation between starbursts and AGN (Taniguchi 2004). However, it is still a matter of debate whether the AGN activity is triggering and causing the starbursts or that the AGN is being fuelled by these massive star forming regions. One thing is certain; there is a growing number of galaxies from different samples that exhibit simultaneous starburst and AGN activity. If there is a causal relation between them, then the question is with which trend. The evolution of supermassive black holes, AGN feeding and feedback to the interstellar medium, as well as the role played by the environment for the formation of stars, are all relevant issues in the physics of starbursts.

### 1.5.3 Feedback: radiative and mechanical

Accreting black holes and massive stars in dense stellar populations produce a large amount of energy. During their active episodes they can emit radiation, produce shock waves, and chemically enrich their surroundings. A lot of this energy is deposited back into the environment from where it was generated. Feedback is a phenomenon where the energy that is created within a system, is fed back into the system itself, altering the ambient conditions and, thereby, influencing the occurrences of the same phenomenon in the present or future. Two types of feedback, radiative and mechanical, are very important in active galaxies during the evolution of star forming clouds. Each type of feedback can have positive or negative influences on star formation.

Radiative feedback will strongly dominate in the form of X-rays, UV, or cosmic rays. Lower energy radiation, like optical or infrared, despite having a strong presence, will be less important in the thermal balance and the chemistry of an interstellar cloud. This is because of their lower energies and their increased attenuation in a dusty and cloudy environment, although this does add a bit to dust heating. Radiative feedback will primarily heat the system as energy is injected into it. However, even then, it is possible to find new ways to enhance cooling. Radiation with high energies $E \geq 1$ keV, like X-rays, will ionize atoms. Ions are much more reactive than neutral atoms and can easily form molecules with other elements. As such, new and more molecules will form, helping the system to cool. On the other hand, radiation with lower energies $E \sim 5 - 100$ eV, i.e., UV, is more destructive to molecules because of large photo-dissociation cross sections.

The strongest X-rays are produced in accretion disks of black holes. A $10^7$ solar mass black hole accreting at 100% Eddington is able to produce a flux of 100 erg $s^{-1}$ $cm^{-2}$ at a distance of 100 pc (Meijerink et al. 2007). X-rays can dominate the thermal balance in AGN upto column densities of $10^{24}$ $cm^{-2}$ and distances of 300 pc (Schleicher et al. 2010b). These regions are known as X-ray dominated regions (XDRs, Lepp & Dalgarno 1996; Maloney et al. 1996), and heating is dominated by photo-ionization. UV radiation has a much smaller penetration depth, which can go upto column densities of $10^{22}$ $cm^{-2}$ (Meijerink & Spaans 2005), at solar metallicity. Therefore, it is most effective if the radiation source is nearby. So UV radiation mainly dominates the chemistry of cloud surfaces. UV radiation heats the gas up to a few thousand K through photo-electric emission from (small) dust grains. The regions where UV radiation dominates, are called photon dominated regions (PDRs, Hollenbach & Tielens 1999). Cosmic rays are not a form of electromagnetic radiation. They are rather energetic charged subatomic particles, mostly protons, moving at relativistic speeds. Cosmic rays can pierce through very large columns of gas and transfer their energy through collisions. Their effect is more subtle and they heat the gas in a molecular cloud almost uniformally. As such, they set the minimum attainable temperature in these systems (Goldsmith & Langer 1978; Bergin & Tafalla 2007).

Mechanical feedback can come from shock waves as well as a strong gravitational potential of a nearby black hole or the deep potential wells in starbursts. Gravitational stresses produce shearing motions that enhance the turbulence of a nearby cloud if it does not tear it apart. This also has major consequences for the accretion rates of proto-stars. Proto-stars in deep potential wells are able to accrete more material than their neighbours (Bonnell et al. 2001; Clark et al. 2008a). A contracting, fragmenting molecular cloud, with massive stars in its vicinity, may enjoy strong shock waves as massive stars go supernova. Shock waves can either compress a cloud and trigger collapse or blow material away.

Nuclei of active galaxies, e.g., ULIRGs like Arp 220 and Markarian 231, enjoy these extreme conditions, see van der Werf et al. (2010). Strong feedback effects thus take place there, but are difficult to observe directly. Little is known on the IMF in active galaxies, so theory and simulations are needed to guide our understanding.

## 1.6 Numerical simulations

Simulation and modelling is an integral part of scientific study. Numerical simulations are a powerful tool that can help to better understand the behaviour of processes. The need for numerical simulations arises when one is limited in performing a study by conventional means, like through analytical work.

In recent years, a large number of studies have been performed on the formation of stars using numerical simulations (Abel et al. 2000; Omukai & Palla 2001; Klessen 2001; Klessen et al. 2005a, 2007; Bonnell & Rice 2008; Wada 2008; Wada et al. 2009; Bate 2010; Hsu et al. 2010; Krumholz et al. 2010; Girichidis et al. 2011; Pérez-Beaupuits et al. 2011; Clark et al. 2011; Latif et al. 2011; Aykutalp & Spaans 2011a). All these studies address a different aspect in the field or improve on an earlier study. The latter can be done through either adding new physics, more resolution and precision, or by implementing a greater dynamical range.

### 1.6.1 Grid codes

Grid based codes are Eulerian codes where one follows the fluid in the lab-frame. The grid is divided into the requested number of cells and the maximum resolution is based on the smallest cell size. The AMR method is a technique which refines the grid only in the regions of interest, according to a predefined refinement criterion, to minimize computational demand while keeping the resolution high. In this way, one can achieve very high resolution at any required location. One of the major drawbacks of AMR codes is its diffusive nature. Especially when advecting over large numbers of grid cells, numerical diffusion is unavoidable. Higher order interpolation schemes reduce the effect, but the issue remains. The only way to minimize this is to increase resolution. On the other hand, a great strength of grid codes is that they can handle shocks and contact discontinuities very well.

Eulerian codes repetitively solve the mass (continuity), the momentum, and the energy equation. In differential form, these equations, coupled with the Poisson equation for gravity, are given by

$$\frac{\partial\rho}{\partial t} + \nabla\cdot(\rho v) = 0 \tag{1.28}$$

$$\frac{\partial\rho v}{\partial t} + \nabla\cdot(\rho v v) + \nabla P = \rho g \tag{1.29}$$

$$\frac{\partial E}{\partial t} + \nabla\cdot(\mathbf{v}(E+P)) = \rho v\cdot g, \tag{1.30}$$

where E is

$$E = \frac{1}{2}\rho v^2 + \rho\epsilon_{\rm int}, \tag{1.31}$$

and the Poisson equation is defined as

$$\nabla^2\phi = 4\pi G\rho \quad \Rightarrow \quad g = -\nabla\phi. \tag{1.32}$$

In these equations, $\rho$ is the fluid (mass) density, v is the fluid velocity, P is the pressure, $\phi$ is the gravitational potential, $\epsilon_{\rm int}$ is the internal energy per unit mass, and g is the gravitational acceleration.

There is a great diversity of simulation codes available within the astronomical community. Ranging from very specific codes created just to serve the purpose of the author to huge, general purpose numerical codes written by groups of people. FLASH, a grid code designed by Fryxell et al. (2000); Dubey et al. (2009), is an example of a large scale, multiphysics simulation code with a wide international user base. The FLASH code forms the basis of the research done in this thesis and is extended with additional physics in each of the chapters.

## 1.7 Thesis outline

The theory of star formation is an interesting, illustrious subject with a lot of discoveries still ahead. In this thesis, the effects of environmental influences on the formation of stars are studied and the results analyzed. The main focus lies on the initial mass function of stars, specifically the formation of stars in extreme environments. Each chapter focuses on a different aspect of star formation and each tells its own tale.

**Chapter 2:**
In the second chapter of this thesis, I focus on the evolution of a giant, r = 10 pc, molecular cloud in a metal-deficient environment. I investigate how metallicity (Z) and an initial rotational moment ($\beta$) affects the fragmentation of this molecular cloud into smaller, but denser, pre-stellar cores. The dependence of molecular cloud fragmentation is tested on the ambient conditions as they pertain to starburst and dwarf galaxy regions, which are then compared against the well known conditions of the

Milky Way. To properly treat the thermal balance, I use a cooling function, created by Meijerink & Spaans (2005), that strongly depends on metallicity. The effects of dust and cosmic rays are also included in the calculations. I simulate the collapsing cloud with these cooling functions for four different metallicities $Z/Z_{\odot} = 1, 10^{-1}$, $10^{-2}$, $10^{-3}$ and for each metallicity condition I consider five rotational energies, i.e., $\beta = 10^{-1}, 10^{-2}, 10^{-3}, 10^{-4}, 0$, where $\beta$ is the initial ratio of rotational to gravitational energy.

**Chapter 3:**
In the third chapter of this thesis, I study a smaller, r = 0.33 pc, turbulent molecular core which exists in the vicinity, at d = 10 pc, of an active black hole. I investigate the effects when this cloud core is irradiated by X-rays, with a flux of 160 erg $s^{-1}$ $cm^{-2}$, emanating from the accretion disk of the black hole. The main question of this exercise is whether star formation is significantly affected by hard X-rays. I perform a full radiative transfer calculation to obtain the proper temperatures inside the molecular cloud by using an XDR code (Meijerink & Spaans 2005). In this study, my focus lies on the emerging IMF in an X-ray dominated region and I assess whether it deviates from a Salpeter shape.

**Chapter 4:**
In the fourth chapter of this thesis, I continue my work on molecular cores in AGN. This chapter directly follows the previous chapter but expands the study on the feedback effects in extreme environments. I now also incorporate the effects of gravitational shear (mechanical feedback), cosmic rays, UV, and varying X-ray fluxes (radiative feedback). A parameter study of 42 different 3D hydrodynamical simulations is performed in order to capture the qualitative and the quantitative effects that the environment imparts on interstellar clouds inside active galaxies. I look at the change in the equation of state and its role in the dynamics of the cloud core. I also analyze how the phase-diagrams, the star formation efficiencies, and the initial mass functions are influenced in active galactic environments.

**Chapter 5:**
In the fifth chapter of this thesis, I combine the results of the previous chapters and evaluate them from a general perspective. I summarize my main findings and give my best answer to the principal question of this work: what is the IMF in active galactic environments? I take the opportunity to discuss what other research paths can be taken to study the origin of, and variations in, the IMF, specifically the role of magnetic fields.